ON THE EVOLUTION OF THE LIGHT ELEMENTS
I. D, ^{3}He, AND ^{4}He

Brian D. Fields

Department of Physics
The University of Chicago, Chicago, IL 60637-1433

and

Department of Physics
The University of Notre Dame
Notre Dame, IN 46556

ABSTRACT

The light elements D, ^{3}He, ^{4}He, and ^{7}Li are produced in big bang nucleosynthesis and undergo changes in their abundances due to galactic processing. Since one may observe most of these elements only in contemporary environments, knowledge of the intervening evolution is necessary for determining the observational constraints on primordial nucleosynthesis. Chemical and stellar evolution model dependences in light element evolution are systematically investigated via a comparison of 1460 possible chemical evolution scenarios and of stellar nucleosynthesis yields, all of which have been selected to fit solar neighborhood C, N, O, and Fe abundances as well as the observed local gas density and gas mass fraction. The light element evolution and solar system yields in these models are found to span a wide range, explicitly demonstrating the model dependence. The range of model dependence for D, ^{3}He, and ^{4}He solar abundances is calculated, and its sensitivity to the heavy element constraints is noted. The chemical evolution contribution to the uncertainty in the observed primordial light element abundances is estimated, and the effects of this uncertainty on big bang nucleosynthesis results are discussed. The predictions for the light elements are found to be correlated; the extent and physical origin of these correlations is discussed. D and ^{3}He evolution is found to have significant model dependence, however, the dominant factor determining their solar and interstellar abundances is their primordial abundance. In addition, ^{3}He is found to be very sensitive to the details of processing in low mass stars. ^{4}He yields are shown to be very model dependent; in particular, both the introduction of mass loss and the possibly very high ^{4}He stellar yields in the poorly understood mass range of $\sim 8 - 12 \ M_\odot$ can lead to large enhancement of ^{4}He production and can lead to large slopes of $\Delta Y/\Delta N$ and $\Delta Y/\Delta O$. It is found that the inclusion of secondary nitrogen leads to only a small distortion in the low metallicity $Y - $ N relation if there is also a significant contribution of primary N, as required by observations.

Subject headings: Galaxies:evolution—nuclear reactions, nucelosynthesis, abundances

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1. Introduction

A central element of the hot big bang cosmological framework is the prediction of light element (D, ^{3}He, ^{4}He, and ^{7}Li) synthesis during the epoch of big bang nucleosynthesis (hereafter BBN). This prediction is tested against observations of the light elements abundances, and the agreement between theory and observation provides major empirical support for the big bang. The comparison between prediction and measurement is not immediate, however. Many of the observations are of environments that are contemporary to the present, or nearly so, and galactic astrophysical production and destruction of the light elements may have changed their abundances. An understanding of these galactic processes is thus needed to in order to accurately interpret the light element observations and so to realistically constrain BBN. The study of the nuclear history of galaxies is the domain of galactic chemical evolution; here we will re-examine the galactic chemical evolution of the light elements.

Models for galactic chemical evolution were first constructed in the 1970's and for the most part studied the nucleosynthesis history of the heavy elements. Since such pioneering work as Cameron & Truran (1971), Talbot & Arnett (1971), and Tinsley (1980, 1972), chemical evolution has come to do a surprisingly good job of reproducing many solar abundance patterns and even many abundance histories as observed in and around the solar neighborhood (see, e.g., the recent work of Timmes, Woosley, & Weaver 1995; Matteucci & Francois 1989; and Ferrini et al. 1992). These models also give reasonable predictions for stellar luminosity evolution (see, e.g. Tinsley 1980; Colin & Schramm 1993), and underpin the study of nucleocosmochronology (e.g. Thielemann et al. 1993).

The chemical evolution of the light elements has been studied since the work of Cameron & Truran (1971), and has been used to determine primordial abundances since the work of Reeves, Audouze, Fowler, & Schramm (1973) and Audouze & Tinsley (1974). These seminal papers made a critical comparison of light element abundances from deuterium to boron with the various cosmological and astrophysical processes suggested for their production. Since this early work there have been many calculations of light element chemical evolution, usually focusing on a few nuclides (see references in §§2.3.–2.5.1.). In this paper we will again take a comprehensive look at all of the light elements to see how consistent a picture of their joint chemical evolution can be made.

Chemical evolution models are physically based on important mass conservation principles, and have been refined considerably over the years. Unfortunately, though, chemical evolution is still a somewhat ambiguous enterprise. One should always be mindful of the uncertainties in the conventional prescription for these calculations. These uncertainties stem from unknown character of star formation rates and mass distributions, galactic inflow and outflow of gas, and from unresolved model dependent effects in stellar evolution nucleosynthesis calculations. Lacking a first principles calculation of these phenomena, we are forced to model them in a phenomenological or schematic way.

As a result of these uncertainties, the commonly adopted framework for chemical evolution contains features that are not well understood in a theoretical and/or empirical way. Thus within this framework, there are a great many ways to construct specific models, and work has been done using a diverse variety of the possible scenarios. Encouragingly, the overall framework has been quite successful despite these difficulties. And indeed chemical evolution is a always a part of comparison of nucleosynthesis calculations with data (though sometimes its role is only implicit).

The effects of chemical evolution model dependences have been investigated to varying degrees. However, the larger studies of these effects

(e.g. that of Tosi 1988) did not concentrate on the light elements, and work on light element evolution has not examined model uncertainties as broadly. There has not yet been an emphasis on the full range of uncertainty in a comprehensive description of light element chemical evolution.

In this paper we construct a chemical evolution framework for the evolution of all of lthe light elements, and we obtain estimates of the range of model uncertainties. We will follow the history of all of the light elements, D through ^{11}B. We also include the non-cosmological "heavy" elements ^{12}C, ^{14}N, ^{16}O, and ^{56}Fe, which serve as tests of potential models. We systematically examine a wide variety of model assumptions, examining many chemical evolution prescriptions and parameterizations (in the spirit of Tosi 1988), as well as including many suggested galactic production and destruction mechanisms for each of the light elements (in the spirit of Reeves, Audouze, Fowler, & Schramm 1973).

Our strategy is to choose a broad chemical evolution framework, allowing for many features that authors have considered in examining the histories of specific light elements. These features are chosen to be inclusive, without regard for their compatibility. With these we form a very large (~ 1200) set of models which include all combinations of these features. We then constrain these models to fit observed solar neighborhood characteristics, apart from those of the light elements. This leaves ~ 120 potential models which are viable descriptions of solar neighborhood chemical evolution (again, light element behavior aside). To each of these successful models we add a range of 18 variants of light element evolution. We then examine the predictions of this large battery of light element evolutionary schemes. Specifically: (1) we see which models fit light element observations; (2) we estimate the chemical evolution uncertainty and discuss its influence on the use of BBN observational constraints; and finally (3) we note correlations between the evolution of different light element

nuclides. As part of this analysis we will consider as well the uncertainties in the solar system constraints themselves.

In addition to these considerations of the whole set of the light elements, we will also address specific issues particular to each of them. For example, recent observations of high-redshift quasar absorption line systems have approached the precision needed to give meaningful D abundances. This is an exciting prospect indeed, as these high-z environments are much more pristine than contemporary ones, and so give a new window to a much earlier epoch in D evolution (though still $\gtrsim 1$ Gyr after BBN). The initial reported abundances by Songalia, Cowie, Hogan, & Rugers (1994) and by Carswell, Rauch, Weymann, Cooke, & Webb (1994) were tantalizingly high if taken at face value (see, e.g. Kernan & Krauss 1994; Cassé & Vangioni-Flam 1994; Steigman 1994). However, the observers themselves have been careful to caution that the absorption line systems may be subject to contamination, and so the derived D abundances are best understood as upper limits to the actual levels of D/H, a caution strengthed by recent reports of a low D abundance in a different line of sight (Tytler & Fan 1994). Clearly, the present situation is in flux and more observations are needed to determine the D abundance in these systems accurately; but just as clearly, this technique is quite promising and holds the key for a very accurate D abundances at very early epochs. We will examine D evolution in detail in this paper, and comment on the role of it chemical evolution in interpreting the quasar absorption line system observational program.

These new constraints on D evolution also call attention to conventional accounts of the evolution of ^{3}He. Recent improved observations of ^{3}He in Galactic H II regions can be puzzling in light of preceding discussions of this element. The data suggest that ^{3}He shows spatial variations, and that the abundance of ^{3}He today in the ISM is not significantly larger than that at the forma-

tion of the solar system—indeed, in some places, ^{3}He is even seen in (possibly) smaller amounts. This may call into question (Hogan 1995) the conventional wisdom about ^{3}He production in low to intermediate mass stars (Iben & Truran 1978). On the other hand, the high ^{3}He abundance seen in planetary nebulae (Rood, Bania, & Wilson 1992) seems to support the notion of production in low mass stars at a level roughly consistent with that predicted by Iben & Truran 1978. Thus the H II region data may more likely point to a need to re-examine ^{3}He evolution in high mass stars (Olive, Rood, Schramm, Truran, & Vangioni-Flam 1994). At any rate, the issue of low mass stellar production of ^{3}He is crucial and we will address it explicitly.

The much more abundant helium isotope, ^{4}He, is important because it is so powerful. Notably, an accurate assay of the ^{4}He content of the universe provides a bound on N_ν, the number of light neutrino families, as first shown by Steigman, Schramm, and Gunn (1977). Thus there are continual improvements in the accuracy of theoretical predictions for ^{4}He (e.g. Dicus et al. 1982; Walker et al. 1991; Kernan 1993; Seckel 1994; Fields, Dodelson, & Turner 1993), as well as in the observed abundances (Skillman et al. 1994). A pressing issue now is the systematic uncertainty in these measurements (e.g. Copi, Schramm, & Turner 1995; Sasselov & Goldwirth 1995). There is also discussion of the correct means of extrapolating the primordial ^{4}He abundance, an issue in which chemical evolution plays a crucial role.

In this paper we will describe our framework for constructing models of light element chemical evolution, and summarize our overall results on model dependences. We will then discuss in more detail our results for D, ^{3}He, and ^{4}He evolution and its uncertainties. In a forthcoming work (Fields 1995, hereafter Paper II) we will discuss our results for Li, Be, and B evolution, with particular focus on the range of available models for cosmic ray nucleosynthesis of these elements.

2. Light Element Evolution

2.1. BBN Predictions

The BBN predictions are crucial to our considerations, as the primordial yields set the initial conditions for the chemical evolution calculation. The basic BBN calculation itself has not changed dramatically since the calculation of Wagoner, Fowler, & Hoyle (1967). For reviews we refer the reader to, e.g., Schramm & Wagoner (1977), Boesgaard & Steigman (1985), and to Smith, Kawano, & Malaney (1993). For recent calculations see Walker et al. (1991), Krauss & Kernan (1994), and Copi, Schramm & Turner (1995).

We consider only the simplest and "standard" case, namely that of BBN in a universe that is spatially homogeneous in all particle species. Some models with baryon inhomogeneities are still allowed by the data, but even these more complicated cases lead to the same basic conclusions, as discussed in, e.g., Malaney & Mathews (1993) and Thomas et al. (1993).

In the standard calculation, the only free parameter is $\eta \equiv n_B/n_\gamma$, the baryon-to-photon ratio. In Figure 1 we plot the light element abundances as a function of η; for reference, Copi, Schramm, & Turner (1995) found concordance between theory and data for the range

$$2.4 \times 10^{-10} \leq \eta \leq 6 \times 10^{-10} \qquad (1)$$

which we have indicated on the plot.

For our purposes there are several salient features to notice. First, the ^{4}He abundance does not depend strongly on η, with variations only at the 10% level; however, we will be interested in the ^{4}He abundance to this level. Even more evident is the decline in D and ^{3}He abundance with increasing η, allowing a large range of initial values, particularly for D which varies more strongly with η. This point is important, as we will find that the a requirement of significant depletion of initially high D down to the solar and interstellar abundance is a serious constraint on chemical evolution models.

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As we will see, the combined abundance D+^{3}He is a most useful constraint on η. Note that for the concordant range of η that this sum is dominated by the contribution from D. Knowledge of the primordial level of D is tantamount to knowledge of D+^{3}He; thus the D abundances obtained from quasar absorption line systems provides a constraint on this sum. Thus an accurate quasar measurement would provide a crucial complement to the solar system and interstellar D and ^{3}He abundances.

2.2. Chemical Evolution Framework

The basic theoretical framework and observational constraints on models of chemical evolution is reviewed in Tinsley (1980). We will, for the most part, adopt her notation. Our goal is to trace the history of the Galaxy's isotopic content. Ideally, we wish to model formation of stars in inhomogeneously distributed gas clouds, their return of material during their lifetime to the ISM via winds, and finally their deaths, which occur in stochastic bursts and return processed and unprocessed material to the ISM. In principle, all of these processes can depend on many variables, e.g. the metallicity of the gas, the temperature of the gas, and the strength and configuration of the galactic magnetic fields, to name a few.

However, this depiction is not computationally feasible. To reduce to problem to a tractable one, we will employ several simplifying assumptions: (1) the Galaxy (or at least the solar neighborhood) is spatially homogeneous; (2) stars are born according to a given mass distribution, and their lifetimes are a function of mass only; and (3) all of the stellar ejecta are returned at once at the time of the star's death, and instantaneously mixed in the interstellar medium (ISM). These assumptions, though strong, are adequate for many applications, and in any case are necessary to proceed.

We will employ the standard convention for the stellar creation rate C, which gives the number of stars born (per pc^2) per Gyr, in dividing

it into a star formation rate (hereafter SFR, ψ) and an initial mass function (hereafter IMF, ϕ)

$$C(m,t)\,dm\,dt = \psi(t)\phi(m)\,dm\,dt \ . \qquad (2)$$

The SFR has units of $[M_\odot\,\mathrm{Gyr}^{-1}\mathrm{pc}^{-1}]$, and gives the mass of ISM material going into stars per unit time. The IMF gives the mass distribution with which these stars are formed. Questions of the time constancy of IMF have been to answer theoretically or observationally (see, e.g. Scalo 1986), although recent work (Beers, Preston, & Shectman 1992) supports the conventional assumption of its constancy, which we will follow.

The literature is fraught with unfortunate variations in the convention for IMF normalization; we follow Tinsley (1980) in choosing

$$\int_{m_{\mathrm{low}}}^{m_{\mathrm{up}}} dm\,m\,\phi(m) \equiv 1 \qquad (3)$$

The lower and upper mass limits, m_{low} and m_{up}, are not well constrained by data, and so will be parameters we will have to choose.

We will use a one zone model for the solar neighborhood, treating it as homogeneous and instantaneously mixed. We will track the total disk mass in terms of its surface mass density σ_{tot}, which evolves by

$$\frac{d\sigma_{\mathrm{tot}}}{dt} = f \qquad (4)$$

where f is the rate for the infall of new material due to, e.g. the collapse of the disk or to accretion of extragalactic material, in units $[M_\odot\,\mathrm{Gyr}^{-1}\,\mathrm{pc}^{-2}]$.

We put the basic constituents of matter as stars (including their remnants) and gas,

$$\sigma_{\mathrm{tot}} = \sigma_{\mathrm{s}} + \sigma_{\mathrm{g}} \ . \qquad (5)$$

Of these, we will explicitly track the gas, which evolves according to

$$\frac{d\sigma_{\mathrm{g}}}{dt} = -\psi + E + f \qquad (6)$$

where

$$E(t) = \int_{m_t}^{m_{\mathrm{up}}} dm\ m^{\mathrm{ej}}\ \phi(m)\ \psi\left(t - \tau_m\right) \qquad (7)$$

is the rate at which dying stars eject processed material back into the ISM. The lower mass limit, m_t, is the stellar mass m for which $\tau_m = t$, i.e. the smallest mass star which dies just at time t after being created at $t = 0$.

We will define the mass fraction

$$X_i \equiv \frac{\sigma_i}{\sigma_{\mathrm{tot}}}\ , \qquad \sum_i X_i \equiv 1 \qquad (8)$$

and we write

$$\frac{d\left(\sigma_{\mathrm{g}} X_i\right)}{dt} = -\psi X_i + E_i + f X_i^{\mathrm{inf}} \qquad (9)$$

where

$$E_i(t) = \int_{m_t}^{m_{\mathrm{up}}} dm\ m_i^{\mathrm{ej}}\phi(m)\psi\left(t - \tau_m\right) \qquad (10)$$

is the rate of the ejection of isotope i to the ISM.

Note that the presence of the "retarded time" $t - \tau_m$ makes the equations unwieldy. In practice one often makes ψ a function of σ_{g} and/or σ_{tot}, thus rendering the basic equations as integro-differential and hence without an analytic solution We have therefore numerically implemented the preceding equations, taking into account their full integro-differential nature. Our results are all based on these calculations.

For comparison with calculations using the instantaneous recycling approximation (IRA), it is useful to define the return fraction

$$R = \int_{m_{\mathrm{low}}}^{m_{\mathrm{up}}} dm\ m^{\mathrm{ej}}\phi(m) \qquad (11)$$

the fraction of mass going into a generation of stars that will eventually be ejected back into the ISM. Note that $R < 1$ and that it is very sensitive to the choice of IMF form and mass limits. We will also use $\mu \equiv \sigma_{\mathrm{g}}/\sigma_{\mathrm{tot}}$ for the gas mass fraction.

Of the model features we have described, those most influential on the abundance predictions are the IMF, the SFR, infall, and the adopted nucleosynthesis yields. Of these, the IMF, SFR, and infall (or outflow) are not well understood theoretically, and so one must select a phenomenological prescription for each. The stellar nucleosynthesis inputs come from model calculations for different stellar masses (and sometimes different metallicities). For the most part only a limited number of metallicities are used, thus forcing chemical evolution models to interpolate and extrapolate. A better treatment would iteratively run both stellar models and chemical evolution models; such a program has recently been done by Timmes, Woosley, & Weaver (1995), who emphasized heavy element yields which they calculated self-consistently using a supernova code.

Of the other input parameters, two are cosmological. The first is of course the BBN result of the initial abundances; our approach to this is discussed below (§4.). The other cosmological parameter is the age of Galaxy including the collapse of the disk; we will take this 15 Gyr, and our results are not very sensitive to this choice. We will take the age of the earth to be 4.6 Gyr. Finally, we will use the τ_m vs. m relation of Scalo (1986).

Having selected a set of models representing the range of choices of the major model features—IMF, SFR, and nucleosynthesis—we will want to constrain this set with observational data. Specifically, we will test these against solar system abundances of CNOFe, as well as the gas mass fraction. The confrontation with observations will select a set of viable models which we will then use to test different models of light element abundances. Thus we will want to include in our initial suite of models all variations of the major features— the IMF, SFR, and nucleosynthesis—which have been proposed to be well-suited to fit individual light elements. Therefore, we now will review the chemical evolution of the light elements, to identify features to include in the initial set of chemical evolution features.

2.3. Deuterium

2.3.1. D Data

Traditionally, D has been measured in the solar system and in the ISM. Important evidence for the solar system D abundance came as a result of the Apollo 11 mission (Geiss & Reeves 1972), giving a measurement of the D+^{3}He content in the solar wind. These data are complemented by measurements of meteorites (Black 1972), which allows determination of the presolar D level. The best numbers now are that the presolar D abundance is $(D/H)_\odot = 2.6 \pm 1.0$ (Geiss 1993).

The D abundance in the ISM was first measured by Rogerson & York (1973). It has most recently been observed by the *Hubble Space Telescope*, with Linsky et al. (1993) reporting $(D/H)_{ISM} = 1.5^{+0.07}_{-0.18}$. This single, remarkably precise measurement will prove to set strong constraints on chemical evolution models.

Most D measurements to date have been of local and relatively recent cosmological epochs. However, exciting new observations have reported possible detection of D in quasar absorption line systems (QSOALS). These systems considered lie at high redshift ($z > 3$), and so are very young (1–3 Gyr). Thus one expects D to have suffered much less depletion. Indeed, initial reports (Songalia, Cowie, Hogan, & Rugers 1994; Carswell, Rauch, Weymann, Cooke, & Webb 1994) suggested the possibility of a very high D, at levels of $D/H \lesssim 2.5 \times 10^{-4}$. Both of these groups have been careful to caution that this result can only be regarded as an upper bound on D, as low column density in intervening clouds can contribute significantly to the putative D line. Indeed, this may be the case for these measurements, as Tytler & Fan (1994) have recently observed a different line of sight and in a high-redshift absorption system claim a preliminary D abundance of $D/H \simeq 2 \times 10^{-5}$. The situation may be more muddled still, as it has recently been suggested (Levshakov & Takahara 1995) that im-

proper modeling of the turbulent characteristics of these systems could lead to very large erros in the derived D abundance. We will thus take the prudent route of not using the few D abundances obtained in this manner, and instead concentrate on better-determined the solar and interstellar values.

2.3.2. D Chemical Evolution

Deuterium was originally thought not to be primordial, but to be produced in the T-Tauri phase of stellar evolution (Fowler, Greenstein, & Hoyle 1962). However, this process was shown to fail (Ryter, Reeves, Gradstajn, & Audouze 1970), and Reeves, Audouze, Fowler, & Schramm (1973) argued that D was likely of cosmological origin. This argument was cemented by the work of Epstein, Lattimer, & Schramm (1976), who showed that not only does stellar burning destroy D, but *any* high energy astrophysical process destroys it as well—except the big bang.

Indeed, as recently emphasized by Copi, Schramm, & Turner (1995), D is best baryometer of BBN in that the constraints on η which arise from D observations are particularly reliable. This reliability stems from the uniquely straightforward evolution of D in galaxies: all of the D in the universe was produced in BBN, and all subsequent processes destroy it. This fundamental fact greatly simplifies the chemical evolution of D, which has no sources and so is absent from any processed material that is returned to the ISM. The D abundance thus provides a clean indication of the amount of material that has not been processed in stars. (The galactic evolution of D ^{3}He, and ^{4}He is summarized in table 1).

To get a feel for D evolution, we note that in the instantaneous recycling approximation, the D evolution can be solved exactly as

$$\frac{X_2}{X_2^p} = \mu^{-R/(1-R)} \quad . \tag{12}$$

Note that this result does not depend on the form of the star formation rate ψ. Furthermore, since

the the gas fraction μ is constrained by observation, the entire character of the depletion comes from R, the return fraction. This is in turn completely determined by the initial mass function. Physically, the D abundance is a measure of the fraction of gas which remains unprocessed. Thus as material is cycled into stars, the D depletion will depend on how much (D-free) material is returned to the ISM to dilute the unprocessed component.

The main chemical evolution feature to consider in order to capture the range of D evolution is that which determines the return fraction R, namely the IMF. We will want to consider the gamut of possible IMF forms and mass limits. Indeed, we will expect all the light element evolution to depend strongly on the adopted IMF.

2.4. Helium–3

2.4.1. ^{3}He Data

The solar system information on both ^{3}He and D come together and are derived from measurements of the solar wind and meteorites (Geiss 1993; Black 1972); the presolar abundance is $(^3\text{He/H})_\odot = 1.5 \pm 0.3$. In the ISM, ^{3}He is observed in Galactic H II regions, and measured via its hyperfine line. Although these measurements are difficult, they apparently show real variations, with $(^3\text{He/H})_{\text{ISM}} \sim (1-5) \times 10^{-5}$ (Balser, et al. 1994). ^{3}He has also been detected by Rood, Bania, & Wilson (1992) in a planetary nebula, and is found at very high levels: $^3\text{He/H} \sim 10^{-3}$.

2.4.2. ^{3}He Chemical Evolution

As noted in §2.4., BBN yields a fairly low ^{3}He abundance, but not much lower than contemporary observations; i.e. BBN theory does not leave a lot of room for the increase of ^{3}He. On the other hand, stellar models firmly show that all initial D in a star is processed to ^{3}He in the pre-main sequence phase when the star is fully convective. Some of this ^{3}He survives to be ejected at the star's death, and indeed in low mass stars are

likely to be ^{3}He sources. Thus while D suffers astration and ^{3}He is produced over time, the sum of the two, D+^{3}He, is more stable; furthermore, D+^{3}He is first dominated by D and then ^{3}He increases until roughly equal measures of each comprise the present ISM abundance.

Yang et al. (1984) get a bound on $(\text{D}+^3\text{He})_\text{p}$ by noting that in all stars, any initial D goes to ^{3}He in the pre-main sequence phase. Some fraction $g_3(m) < 1$ of that ^{3}He survives in the ejecta at the star's death. If we consider one stellar generation, and let $\langle g_3 \rangle$ be the average of $g_3(m)$ over the mass function, it may be shown that

$$\left(\frac{\text{D} + {}^3\text{He}}{\text{H}}\right)_\text{p} \leq \frac{1}{\langle g_3 \rangle}\left(\frac{{}^3\text{He}}{\text{H}}\right)_1 + \left(\frac{\text{D}}{\text{H}}\right)_1 \quad (13)$$

(Yang et al 1984). The inequality (eq. 13) comes from disregarding ^{3}He production, and so guaranteeing that X_{23} decreases with time.

We may get another perspective on this point by seeing how the D+^{3}He argument plays out not for a single stellar generation, but in the approximation of instantaneous recycling (see preceding section) for a closed box chemical evolution model. We have

$$\frac{d\left(\sigma_\text{g} X_3\right)}{dt} = -\psi X_3 + R\langle g_3 \rangle \psi \left(\frac{3}{2} X_\text{D} + X_3\right) \quad (14)$$

and when substituting from eq. 9 we have

$$\frac{(\frac{3}{2} X_2 + X_3)}{(\frac{3}{2} X_2 + X_3)_\text{p}} = \left(\frac{X_2}{X_2^\text{p}}\right)^{1-\langle g_3 \rangle} \quad (15)$$

and so in term of observable number ratios,

$$\begin{aligned}
\frac{(\text{D} + {}^3\text{He})}{(\text{D} + {}^3\text{He})_\text{p}} &= \left(\frac{\text{D}}{\text{D}_\text{p}}\right)^{1-\langle g_3 \rangle} \left(\frac{X_1}{X_\text{p}}\right)^{\langle g_3 \rangle} \\
&\approx \left(\frac{\text{D}}{\text{D}_\text{p}}\right)^{1-\langle g_3 \rangle} \quad (16)
\end{aligned}$$

D+^{3}He increases (decreases) with time if $\langle g_3 \rangle > 1$ ($\langle g_3 \rangle < 1$). Note, however, that because low

mass stars are crucial here, that the IRA is inappropriate. At any rate, there is a close relationship between D and ^{3}He, and we see that it is crucial to know whether ^{3}He is created or destroyed with time.

The physical explanation for the different fate of ^{3}He in high and low mass stars is that for $m < 2M_\odot$, $p - p$ is the dominant form of hydrogen burning, and the $p(p, \nu e^+)D(p, \gamma)^3$He chain makes ^{3}He. However, for $m > 2M_\odot$, the CNO cycle dominates the hydrogen burning, and ^{3}He is destroyed. Indeed, Iben & Truran's (1978) calculation for low mass stars suggests that ^{3}He is produced copiously. Dearborn, Schramm, & Steigman (1986) find that ^{3}He is destroyed in high mass stars, particularly at low metallicity. Woosley & Weaver (1994) give more detailed but qualitatively similar results.

These conventional notions about ^{3}He evolution have recently been called into question as the interstellar ^{3}He abundances have become more accurate. Until recently, the production of ^{3}He in low mass stars was not considered in chemical evolution studies, e.g. that of Steigman & Tosi (1992). Vangioni-Flam, Olive, and Prantzos (1994) let the contribution to $\langle g_3 \rangle$ from low mass stars be a free parameter, and used the Dearborn et al. (1986) high mass yields. They found that to fit the data require net destruction by low mass stars, a mechanism for which is suggested by Hogan (1995). Indeed, with the Iben & Truran (1978) and Dearborn et al. (1986) results, Galactic ^{3}He which starts with the espoused BBN value is apparently overproduced in many chemical evolution models (Olive et al 1994; Galli et al. 1994). However, in the face of this apparent need for ^{3}He destruction is the observed prescence of high ^{3}He in planetary nebulae (Rood, Bania, & Wilson 1992) at a level in rough agreement with Iben & Truran's calculation.

Another ^{3}He puzzle is the observed and apparently real dispersion in ^{3}He abundances. Balser et al. (1994) present this dispersion as a ^{3}He abundance gradient, with ^{3}He decreasing towards the galactic center. This trend runs counter to intuition, if ^{3}He is to be formed by low mass stars which are more prevalent towards the galactic center. Moreover, for stars having lifetimes this long, it is hard to see how this gradient could happen at all, as such stars will have traveled a significant fraction of disk in their lifetimes and so should be well spread throughout the Galaxy.

A possible explanation for "what is wrong with ^{3}He" was offered by Olive et al. (1994). They point out that the correlation of ^{3}He with galactocentric distance might be better understood as a correlation with the mass of the H II region. Then the ^{3}He in the higher mass regions will contain more ejecta from high mass stars. In this scenario, one expects ^{3}He to be lower wherever there is more ejecta from massive stars, i.e. more stellar processing. Thus the ^{3}He abundance should go down towards the center of the galaxy, where the density and so the star formation rates are higher. Note that this could also answer the puzzle of the existence of the gradient despite the long lifetime of low mass stars. The point is that there is not a gradient in ^{3}He production (by low mass stars), but instead there is a gradient in destruction by short–lived, high mass stars.

For our chemical evolution models, we will want to use our apparatus to address the important issue of the effect of low mass processing on ^{3}He evolution, allowing for different possibilities as done by Vangioni-Flam, Olive, & Prantzos (1994). We will also include infall, as was done in Steigman & Tosi (1992).

2.5. Helium–4

2.5.1. ^4He Data

Observations of ^{4}He are well reviewed elsewhere, e.g. Pagel (1993), Pagel et al. (1992), Walker et al. (1991), Davidson & Kinman (1985), Skillman & Kennicutt (1993); Skillman et al. (1994); and Copi et al. (1995). Briefly, accurate determinations of ^{4}He are hard to come by. The best observations are of H II regions, where

He should be ionized and accurate abundances might result. Although there have been observations of ^{4}He in Galactic H II regions, systems with lower metallicity are expected to contain less contamination from stellar ejecta. The sites of choice have proven to be metal poor extragalactic H II regions.[1]

The solar system abundance of ^{4}He is $Y_\odot = 0.274 \pm 0.016$ (Anders & Grevesse 1989). ISM abundance determinations are few, and furthermore, the ionization structure of gas regions causes ^{4}He to be systematically underestimated. (Wilson & Rood 1994). Thus, we have no information on ^{4}He in the ISM that is accurate enough to be an important constraint on chemical evolution models.

2.5.2. ^{4}He Chemical Evolution

As the second most abundant nuclide in the universe, next to hydrogen, ^{4}He comprises about a quarter of the universe's baryonic mass. It is also the only nuclide made in significant amounts in stars of a broad range of masses, with the dominant production in stars of middle masses (Renzini & Voli 1981; Iben & Truran 1978). Despite this nearly ubiquitous stellar synthesis, however, stars only make a small ($< 10\%$) contribution to the total helium abundance (Hoyle and Tayler 1964): most ^{4}He is primordial. Nevertheless, to fully test BBN and moreover to use its power to constrain particle physics and/or test conditions in the early universe, one needs to know the abundance of ^{4}He exceedingly well.

Even in low metallicity H II regions, stellar pollution exists; the question is how best to determine it and so deduce the primordial He abundance. To this end Peimbert & Torres-Peimbert (1974, 1976) noted that He production should be correlated with heavier element production. In

H II regions one observes both He and CNO. Following Peimbert & Torres-Peimbert, one deduces the primordial helium mass fraction Y_p from extrapolating the low metallicity end of the Y vs. Z plot, exploiting the relation

$$Y = Y_p + \frac{dY}{dZ}\,Z \qquad (17)$$

Equation (17) should be valid for sufficiently small Z (and dY/dZ a constant function of Z), for which this procedure should succeed in extrapolating Y_p. The challenge for chemical evolution is to determine which Z are sufficiently small, and which "metals" C, N, and/or O are the best surrogates for Z. The challenge for observers is to make precise enough measurements to allow for a meaningful extrapolation.[2]

Analysis using this technique has for the most part avoided the explicit use of detailed chemical evolution calculations. Instead, the approach has been an empirical one: Y and Z are measured for different regions; the Y vs. Z relation is plotted, fit, and extrapolated to get $Y_p = Y(Z = 0)$. Two complications to this procedure immediately arise. First, one does not directly measure the metal fraction Z, and so one uses as a surrogate either oxygen or nitrogen (and carbon—see Steigman, Gallagher, & Schramm 1989). Second, and moreover, one must adopt an expected Y vs. O (and N or C) relation to use in fitting the data and making the extrapolation.

While everyone recognizes that the fits need not be linear, this is the simplest two parameter fit, and so this has been the first to be tried. A glance at the data (see figure 7) suggests that this might not be a bad first guess. Indeed, the Y vs. O/H relation is well described by a linear fit. In contrast, the case of nitrogen as a metallicity

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[1]Jakobsen et al. (1994) have recently reported detection of ^{4}He in a quasar absorption line system. The uncertainties in the measurement render it a qualitative confirmation of BBN, but with more accuracy this type of observation could be extremely powerful.

[2]The difficulties of measuring He and metal abundances to the desired accuracy (Y_p good to the third decimal place!) cannot be overstated. For attempts to correct for some of these errors by culling points from the data set see Pagel (1993), Pagel & Kazlauskas (1992), Pagel, Simonson, Terlevich, & Edmunds, (1992), and Olive & Steigman (1995).

tracer has generated vigorous debate as to the appropriate fitting function. At issue is whether to choose a linear fit, in which the slope $\Delta Y/\Delta N$ is constant, or a nonlinear one, in which $N \propto Y^2$, and thus the slope is not constant and has the possibility of being quite steep for small N, i.e. at early times (Fuller, Boyd, & Kalen 1991; but see Olive, Steigman, & Walker 1991).

Put differently, the question regarding the $Y - N$ relation is, what is the $N - O$ relation? The nonlinear fit for Y vs. N arises by assuming that N production is proportional to the C and O abundance, leading to the "secondary" relationship $N \propto O^2$. We wish to know the relative size of the primary and secondary stellar contributions to N, an issue not yet resolved. Consequently, we will want to test effects of both primary and secondary N.

Most ^{4}He observations are determined from extragalactic H II regions. To completely understand these abundances would require separate evolutionary models for each observed galaxy. However, we will be fitting our model to the solar system, and so we should not necessarily expect a close fit to the extragalactic data. However, the few models that do exist, e.g. those of Mathews Boyd & Fuller (1993), and Balbes, Boyd, & Mathews (1993), are qualitatively similar to those studied here.

Another outstanding problem in ^{4}He evolution is the theoretical reproduction of the slope $\Delta Y/\Delta Z$. The theoretical dY/dZ depends on how many high mass stars are included in the calculation (e.g. Maeder 1983, 1992, 1993; Brown & Bethe 1994; and Prantzos 1994). Stellar yields for massive stars give a much lower ratio $dY/dZ \sim 1.3$ than that for intermediate and low mass stars, $dY/dZ \sim 6$. Averaged over a typical IMF with mass limits of 0.4 and 100 $M_\odot$ gives

$$dY/dZ \sim 2 \qquad (18)$$

at most (Maeder 1992, 1993). This theoretical slope is to be compared with an observed slope (calculated assuming $Z \propto O$) of $dY/dZ \sim 4 - 6$

(Pagel 1993).

One attempt to address this issue notes that many models for high mass stars (e.g. Weaver & Woosley 1993) do not include mass loss. For stars above $\sim 30~M_\odot$ this can have a significant impact on all abundances but particularly ^{4}He, as suggested by the results of Maeder (1992). Indeed, stars above $\sim 30~M_\odot$ are thought to rapidly lose mass until reaching $30M_\odot$; thus one need only compute stellar yields up to this mass. For our purposes, we will explicitly include the possibility of mass loss in high mass stars.

Finally, an unfortunate aspect of stellar evolution is that there is a great deal of uncertainty as to the gross behavior–let alone the nucleosynthesis yields–of stars in the 8–10 $M_\odot$ mass range (further discussed below, §3.1.1.). It has been suggested (Woosley & Weaver 1986) that these stars might dominantly produce helium. This could have a large effect on the net ^{4}He production and so it is worth trying $8 - 10M_\odot$ yields with very large yields helium, to see the sensitivity to this range.

3. The Suite of Models

Our goal is to examine a broad selection of physically plausible scenarios, including at least the variations that have been commonly used in the literature. This approach is inspired by that of Tosi (1988), although her study made infall a central feature whereas we will emphasize other features as well.

We classify model features as "global" in scope for particular to the light elements. We will designate "global" model features to be: (1) those that affect all elements (e.g. choices of SFR or IMF) and the gas consumption; or (2) those that affect the yields of C, N, O, and Fe (CNOFe), i.e. the nuclides we will be following which are not light elements but serve as chronometers or tracers for the light elements. Upon identifying the set of global model variations, we combine all of these features independently to create a

large set of potential models for the solar neighborhood. By systematically investigating every possible combination of our chosen global features, we will be assured to properly estimate the full model variation inherent within the span of model options we have allowed.

We demand that our results conform to the features of the solar neighborhood. By selecting those models which reproduce the solar characteristics, we will develop a suite of acceptable global models. To each of these global models is then added a wide variety of light element evolutionary prescriptions. We then calculate the chemical evolution for every light element variant, obtaining an inclusive measure of the model dependence of light element evolution.

3.1. Selection of Global Model Features

In her study, Tosi (1988) found that all aspects of chemical evolution models affect the absolute abundances. However, she found that the choice of IMF, and of course the selection of nucleosynthesis yields, had the most profound influence, affecting the abundance ratios as well as the absolute abundances. Therefore we will want to be particularly thorough in our consideration of these model features. Our global model properties are summarized in table 2.

3.1.1. Nucleosynthesis Yields

The tabulated nucleosynthesis yields are those of low to intermediate mass stars, with masses $\sim 0.8 M_\odot$ to $\sim 8 M_\odot$, and high mass stars with masses $M > 12 M_\odot$. The former are understood to be stars whose lives progress to the point of planetary nebula ejection and becoming carbon-oxygen white dwarfs; the latter, high mass stars end their lives explosively as type II supernovae. Standard results for this study are those of Renzini & Voli (1981) for intermediate mass stars, and Weaver & Woosley (1993) for high mass stars.

The uncertainties in the yields pointed out in these references. The Renzini & Voli (1981) results, for example, were found to depend on the adopted parameterization of convection in the form of the mixing length. However, this level of detail is not a major source of uncertainty for our purposes. It would be desirable to compare results for different stellar models and yields; unfortunately this is not possible for the low mass stars, as Renzini & Voli (1981) are the only current models with detailed results reported. In addition, for the massive stars, several groups have published massive star yields; however these models have very different degrees of emphasis on their nuclear yields, with Woosley & Weaver (1993) presenting particularly meticulous results.

As motivated in §2.5., we allow for a wide range of N evolution, including both primary and secondary sources for N. We thus will follow the approach of Mathews, Boyd, & Fuller (1993) in writing the mass fraction of ejected N, X^{ej}, as

$$X^{\mathrm{ej}} = X_0^{\mathrm{ej}} \left(\alpha + \beta \frac{X_C(t)}{X_C^\odot} \right) \qquad (19)$$

where X_0 is the usual N yield, $X_C(t)$ is the calculated C mass fraction and $X_C^\odot$ is the solar carbon mass fraction from Anders & Grevesse (1989). The constants α and β are chosen as follows:

$$(\alpha, \beta) = (1, 0) \; ; \; (0.5, 1) \; ; \; (0, 1) \; ; \; (0, 2) \qquad (20)$$

where the first option is the standard case of purely primary N, the second is a mixed case of some primary and some secondary N, and the final two cases are variation on purely secondary N. Note that our procedure differs somewhat from that of Mathews, Boyd, & Fuller (1993) who used O as the seed nucleus for secondary N, instead of C. Carbon is the more appropriate choice (Audouze, Lequeux, & Vigroux 1975; Vigroux, Audouze, & Lequeux 1976; and Dearborn, Tinsley, & Schramm 1978), and we have used it.

3.1.2. Initial Mass Function

The selection of an initial mass function is a crucial feature of any chemical evolution model.

Unfortunately, there is no convincing physical theory of star formation, and so the nature of the IMF and the star formation rate over the history of the Galaxy are poorly understood. It is unclear, for example, whether the IMF has changed with time, what its upper and lower mass limits are (and have been), and whether it depends upon the composition of the gas which will become stars. Furthermore, it is very difficult to untangle the behavior of the IMF and the SFR, which could in fact be inseparable and better left in the form of a creation function, as in eq. (2).

Lacking any good theoretical or observational guidance as to these questions, we will adopt eq. (2), i.e. we will assume that the IMF is constant in time. We will, however, allow for it to take different forms. We will try the classic Salpeter (1955) function,

$$\phi(m) \propto m^{-(1+x)} \qquad (21)$$

and we will allow the index ("slope") x to vary, trying $x = 1, 1.35$, and 1.7. We will also try an IMF ϕ derived from the observed, present-day mass function (PDMF, ϕ^{pd}), as given by, e.g. Scalo (1986). The relationship between the two depends on whether the stellar lifetime τ_m is short enough that some of the stars of mass m have died; specifically,

$$\phi(m) = \begin{cases} \phi^{\mathrm{pd}}(m)\,, & \tau_m \geq t_0 \\ \phi^{\mathrm{pd}}(m)/\int_{t_0-\tau_m}^{t_0} dt\, b(t)\,, & \tau_m < t_0 \end{cases} \qquad (22)$$

where the age of the Galaxy is t_0 and $b(t) = \psi(t)/\langle\psi\rangle$. We remind the reader that this procedure is only practicable when the star formation rate ψ is a given, explicit function of time,[3] because of the role of ψ in eq. (22).

We also will try different mass limits to the IMF. Here we are guided by common choices that

have appeared in the literature. We will wish to have a broad, medium, and narrow range in mass. We choose limis of

$$(m_l, m_u) = (0.2, 100); (0.1, 60); (0.4, 30) \qquad (23)$$

which we feel to be representative.

3.1.3. Star Formation Rate

As the physics of star formation is ill understood, the form of the SFR is not well set by theory, although there have been some attempts to do so. As with the IMF, there are many variants, but we will select what we feel to be typical choices:

$$\psi = \nu\sigma_{\mathrm{tot}}\left(\frac{\sigma_{\mathrm{g}}}{\sigma_{\mathrm{tot}}}\right)^0 = \nu\sigma_{\mathrm{tot}} \qquad (24)$$

$$\psi = \nu\sigma_{\mathrm{tot}}\left(\frac{\sigma_{\mathrm{g}}}{\sigma_{\mathrm{tot}}}\right)^1 = \nu\sigma_{\mathrm{g}} \qquad (25)$$

$$\psi = \nu\sigma_{\mathrm{tot}}\left(\frac{\sigma_{\mathrm{g}}}{\sigma_{\mathrm{tot}}}\right)^2 \qquad (26)$$

$$\psi = a\,\exp(-t/\tau) \qquad (27)$$

where we pick the $\tau = 7.5$, 15 Gyr as the timescale for the exponential case. Note that these always decrease for a closed box ($\sigma_{\mathrm{tot}} = $ const), but that for infall models with accreting gas, these will show a rise at early times.

3.1.4. Infall

There is very little known about the existence, nature, and composition of infall and/or outflow from our Galaxy. Evidence for infall is perhaps given by the observation of the High and Very High Velocity clouds. However, there is certainly a need for infalling matter at early times, if one views infall to disk as just the outflow from the halo as it collapsed. In this case, one imagines a metal-poor (i.e. BBN composition) infall on a short (~ 2 Gyr) timescale. Another justification for infall is more pragmatic: it provides a possible solution to the G-dwarf problem (discussed in §3.2.).

We will adopt models with and without infall, which we parameterize according to the widely

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[3]In fact, other SFRs can accommodate the Scalo IMF through an iterative process (Mathews, Bazan, & Cowan 1992); such a procedure would however be very time consuming to use with our many models, and is not considered here.

used form $f(t) = f_0 \exp(-t/\tau_{\rm inf})$. We will vary the infall strength f_0, and the timescale $\tau_{\rm inf}$. Namely, we will try $\tau_{\rm inf} = 2, 4$ Gyr, and we will use different f_0 such that infall contributes 50% and 99.9% of the total disk mass today. We will also consider "closed box" models with no infall, i.e. for which $f_0 = 0$.

3.1.5. Supernova Rates

We will assume that all supernovae arise either from the core collapse of massive stars (type II) or from accretion onto a CO white dwarf in a binary system (type Ia). All stars above $\sim 12 M_\odot$ are thought to become type II supernovae, whereas the CO progenitors of type Ia supernovae are arise from binary systems of intermediate $(3 - 8\ M_\odot)$ mass. The rates for these events are given by (Matteucci & Greggio 1986), and are of the form

$$R_{\rm Ia}(t) = \lambda \, \mathcal{F}[\phi, \psi, f] \qquad (28)$$

where $\mathcal{F}$ is a functional of the IMF and SFR, as well as f, the primary-to-secondary mass distribution in binaries (and perhaps also the initial separation; e.g., Mathews et al. 1993).

The amplitude of the type Ia rate is controlled by the dimensionless parameter λ which is the probability for binary systems in the appropriate mass range to undergo supernova events. This number is not well understood theoretically and in practice is adjusted to reproduce the present ratio of type II to type Ia events ~ 1; this ratio is proportional to λ but is much larger due to the larger number of low-mass, potential type Ia progenitors. We will examine the sensitivity of our results to this parameter, by choosing two values for λ: a high value of $\lambda = 0.05$ near the value recommended by Matteucci & Greggio (1986), and a low value $\lambda = 0.007$ used by Timmes, Woosley, & Weaver (1995).

The type II nucleosynthesis yields are those discussed in §3.1.1., and are directly included in eq. (9). Type Ia yields of nuclide i are added by putting

$$E_i^{\rm Ia} = M_i^{\rm ej} R_{\rm Ia} \qquad (29)$$

with $M_i^{\rm ej}$ the mass ejected in i in a type Ia event. We use the yields for the W7 model of Thielemann, Nomoto, & Yokoi (1986).

3.2. Constraints on the Global Models

We first constrain the global models to fit the solar neighborhood properties within fairly generous tolerances, excluding for the moment any consideration of the acceptability of the global models' light element yields. The available constraints on the gross features of the models are $\sigma_{\rm g}^1$, μ_1, and the CNOFe abundances.

Note that we want to be not too restrictive on our models, as the data, even the solar system abundances, are likely to be subject to systematic errors. For example, Olive & Schramm (1981) suggest that the solar system may have been formed from a young OB association; as a result its elemental and isotopic content would be a biased indicator of the larger solar neighborhood. Indeed, it seems unwise to dwell on one constraint at the exclusion of the rest. Consequently, we wish to be generous in choosing models with which to check chemical evolution uncertainties in the light elements.

We will demand that all models fit the local gas mass fraction, which is determined to be $\mu = 0.1 - 0.2$ (Rana 1991). Specifically we require $\mu = 0.13 - 0.17$. This fixes the overall mass consumption. Given that the a central feature explicitly built into chemical evolution models is the mass consumption and the conservation of overall mass, this is a most important constraint. We therefore have tuned all of the SFR normalizations to satisfy this constraint.

As for the quantitative means of evaluating a model, the χ^2 analysis of Tosi (1988) is interesting but perhaps premature given that the data is likely to be fraught with systematic errors. Instead, we will make the generous (and implementationally simple) demand that the solar CNOFe

abundances within a ± 0.4 dex[4] range. This is to be compared with the factor of 2 that Timmes, Woosley, & Weaver (1995) allowed themselves in a model tuned to give good results.

We will also constrain some of our models with observed G-dwarf distribution. G-dwarf stars are long lived, with ages comparable to the Galactic age now. Consequently, their provide a record of the integrated star formation history of the galaxy; the metallicity distribution of these stars is as seen in Figure 9. It has long been known that the simplest models of galactic chemical evolution are unable to reproduce the G-dwarf distribution. In particular, such models cannot avoid overproduction of low metallicity stars, whereas the data shows a dearth of G-dwarfs below a metallicity of [Fe/H].

One of the solutions to this classic "G-dwarf problem" is the introduction of infall. In models with infall, low metallicity stars are still formed, but as material is added subsequently, the fraction of low metallicity G-dwarfs becomes acceptably small. We will thus apply a G-dwarf constraint to our models, but only on models with infall. For these we will demand a solution to the most emphasized part of the G-dwarf problem, namely the need for a small number of low metallicity G-dwarfs. We will ask that our models not exceed the observed fraction of such stars at [Fe/H] = -0.8 by more than a factor of 0.4 dex $\simeq 2.5$; this will prove to be a restrictive condition. We will not make the stronger demand that the models fit the full G-dwarf distribution to a high accuracy, as this is not done even in models which allow themselves tuning which we lack.

We will not impose the G-dwarf constraint on models without infall, as it has long been thought (e.g. Tinsley 1980; Pagel 1989) that such closed box models are unable to meet this constraint

without further modification. On the other hand, Mathews & Bazan (1990) show that the G-dwarf problem is somewhat mollified by the inclusion of metallicity dependent stellar ages, and Sommer-Larsen (1991) has noted a selection bias in favor of high metallicity dwarfs. At any rate, we keep the closed box models since their ubiquitous presence in the literature demands that one examine their results, and again, we want to err on the side of inclusiveness in our selection of viable models.

3.3. Light Element Model Features

Having assembled a set of possible models and determined constraints on their gross characteristics, we now turn to the model features we will use to encompass the possibilities for light element evolution. For the most part, these introduce small enough changes in the CNOFe abundances that each model's CNOFe prediction will remain valid despite the modification of light element evolutionary features. An exception to this rule comes from mass loss.

The evolution of all light elements is affected by their initial BBN abundance as set by the choice of η. We use three different values of η, guided by the analysis of Copi, Schramm, & Turner 1995 (cf. eq. 1). We choose central, high, and low values $\eta = (2, 4, 6) \times 10^{-10}$; these values bracket the Copi, Schramm, & Turner's "sensible" limit. The low limit is intentionally chosen to be slightly below their recommended value (but above their "extreme" lower limit), in order to allow comment on recent speculation of the possibility of a low η.

In addition to systematically varying η, we will include the following features for each element:

3.3.1. D

As we have discussed, D evolution is uniquely simple because there are no stellar sources of D; the range in D results thus stems only from the choice of η and from global model features.

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[4]dex = base 10 logarithmic units; i.e. a logarithmic variation by ± 0.4 dex is a linear variation by a factor of $10^{0.4} = 2.5$

3.3.2. ^3He

For high mass ^{3}He yields we used the results of Weaver & Woosley (1994), which span a grid of masses from $\sim 11 - 40 \ M_\odot$ and very complete range of metallicity from $Z = 0$ to $Z = Z_\odot$. The metallicity dependence of these models is strong and thus a crucial feature to include; these results update the work of Dearborn, Schramm, & Steigman (1986).

As noted in §2.4., ^{3}He evolution hinges on the effect of low mass stars on ^{3}He. To investigate this effect we have chosen an approach similar to that of Vangioni-Flam, Olive, & Prantzos (1994), as well as Olive et al. (1994). Namely, we specify the ^{3}He survival fraction g_3 at $m = 1$, 2, and 3 $M_\odot$. We choose to allow for: (1) modest ^{3}He destruction (Vangioni-Flam, Olive, & Prantzos found they must); (2) "break even," with all ^{3}He surviving but no net production; and (3) modest ^{3}He production. Specifically, the three ranges we choose are

$$g_3 \left(1M_\odot, 2M_\odot, 3M_\odot\right) = \begin{cases} (1, 0.7, 0.7) \\ (1, 1, 1) \\ (1.33, 1.33, 1.33) \end{cases} \quad (30)$$

We chose the production to be small (i.e. somewhat less that the Iben & Truran (1978) yields, corrected to include initial D, would suggest) in order to be conservative. It is clear, as pointed out in Olive et al. (1994) that the yields at face value do not work, assuming the conventional ^{3}He high mass yields. In an effort to allow for uncertainties in the Iben & Truran (1978) calculation, we wish to see if even a modest production can work.

3.3.3. ^4He

As discussed in §2.5., calculations of ^{4}He production must allow for its production by stars having a wide range of masses; because of this broad range, ^{4}He is particularly sensitive to systematic problems in nucleosynthesis yields over the whole mass range. One such problems is that tabulations of stellar nucleosynthesis yields usually do not include results for stars in the range of roughly $8 - 12 \ M_\odot$. There is not a firm understanding of the basic nature of these stars' final fate; whether they become white dwarfs or explode as supernovae is a delicate and unresolved question. The nucleosynthesis results for these stars are therefore quite uncertain, and indeed possibly quite different from those of stars either more or less massive. This is a regrettable state of affairs, as a typical IMF places a good deal of weight on the yields from stars in this range. Thus whatever assumptions one adopts about these yields can prove important for chemical evolution, and so we will examine two possible scenarios.

Traditionally, one obtains yields in this mass range by interpolating the results from the closest tabulated masses. Such an interpolation can be viewed as hedging one's bets about the accuracy of estimates for the onset of core collapse, and we will include this case; however, it has also been suggested that stars in this range produce ^{4}He almost exclusively (Woosley & Weaver 1986). We therefore mock up this behavior by assuming that *all* of the ejecta from these stars is in this form.

Other important model uncertainties regarding ^{4}He production arise due to the omission of mass loss in high mass stellar models Consequently, we use the stated Woosley & Weaver (1993) high mass yields, which do not include the effects of mass loss, but we also characterize the effects of mass loss by adding Maeder's (1992) calculation of these. Note that while Maeder also calculated supernova yields which included mass loss, the models were far less detailed and thorough than those of Woosley & Weaver; thus we choose to import only the mass loss addition.

4. Results

4.1. Effect of the Chemical Evolution Model Features

The effect of our global model features are for the most part well documented in the literature; we will discuss only the highlights. For a more complete discussion we refer the reader to, e.g., Talbot & Arnett (1971), Tinsley (1980), and Tosi (1988). Note that unless otherwise indicated, all of the figures in this section are for models which vary only the parameters discussed while keeping the others constant.

A sample of SFR behaviors appears in Figure 2; Figure 2 (a) displays the case of a closed box model. We see that for SFRs of the form $\psi = \nu\sigma_{tot}\mu^n \propto \sigma_g{}^n$, there are progressively higher initial bursts and more rapid declines for progressively higher n. The presence of infall changes the behavior of the SFRs whose form depends upon the total and/or gas masses. As seen in Figure 2 (b), for $\psi = \nu\sigma_{tot}$ the SFR is monotonically increasing with time as the disk mass grows through infall, in contrast to the decaying SFR forms in a closed box. For $\psi = \nu\sigma_g$ (Figure 2 (c)), in the case of infall the SFR initially rises to a peak, but then falls off as the gas is converted to stars and remnants.

The total and gas mass evolution is illustrated in Figure 3. Our SFR boundary conditions and infall prescription fix the σ_{tot} plot as well as the endpoints of the σ_g plot. Note that the spread in σ_g at $t_1 = 15$ Gyr is an indication of this spread in μ_1, which arises due to slight differences in the SFR tuning for different models. The effect of infall on the gas and total mass is clear from Figure 3. Without infall, the total mass is of course constant, and the gas mass monotonically decreases. With infall, the total mass rises rapidly, with the gas mass initially following, then reaching a peak and turning over.

The relation between iron abundance and time (the "age-metallicity relation") is an important one: iron is readily observable whereas t is not, and the maximal binding energy of Fe ensures that its abundance will increase monotonically with time. Thus the iron abundance is traditionally used as a chronometer (although its origin in both type Ia and type II supernovae makes it a somewhat problematic one, with O perhaps being a more suitable choice. See Wheeler, Sneden, & Truran 1989).

In Figure 4 we plot [Fe/H] versus t for all of our SFR options, with data from Edvardsson et al. (1993). In general, we see that the Fe rises very quickly, with most of the production happening at very early times and very little new Fe being added today. In comparing the curves in Figure 4 to the data it is immediately clear that all of the models are viable. There is also a great deal of scatter in the data, indeed, as much as the overall trend. Edvarsson et al. (1993) emphasize that this scatter is real, and is not well explained in most chemical evolution models which invoke, as we have, the assumption of instantaneous mixing of stellar ejecta at a given epoch. We can only expect to reproduce the average trends in the observations, and one should recall that individual stars can show considerable excursions from the average trend.

The evolution of deuterium provides a good example of the effect of different IMFs; the D history for three different IMFs is displayed in Figure 5 (a). The three choices have different return fractions, with $R = 0.32, 0.42$, and 0.57. As expected (cf. eq. 12), the D evolution is very sensitive to this parameter, with the overall depletion at the solar system (i.e. at $t_\odot = 10.4$ Gyr) varying from a factor of about 1.4 to a factor of 2.3. As we will see, this spread accounts for a large portion of the model uncertainty in the D evolution (for a fixed η). Note that the largest D depletion comes from using the Scalo (1986) IMF, which is derived from the present-day mass function. This IMF is bimodal and so has a lot of power in the heaviest stars which eject most of their mass; consequently, this yields a large return fraction and considerable D depletion.

In Figure 5 (b) we show ^{3}He and D+^{3}He for

the same models. The tradeoff between D and ^{3}He is clearly seen, with ^{3}He being progressively higher in models having progressively lower D. Indeed, despite the large variation in the individual D and ^{3}He cases for the three IMF choices, we see that all show a very similar—and nearly constant—evolution of D+^{3}He. In Figure 5 (c) we show the ^{4}He evolution for the same models. These models show a correlation between the D and ^{4}He evolution, as the models with larger D depletion also have larger ^{4}He production. We will investigate this relationship in detail in §4.3..

Deuterium together with ^{3}He provides a good barometer of the effects of infall on the evolution of an isotope. The infall of primordial material enriches the depleted Galactic D, while it dilutes the increasing Galactic ^{3}He. This behavior is evident in Figure 6 which shows the slower decline of D in the infall models, mirrored by the inhibited growth of ^{3}He. Note, however, that in the case of D the spread (at the time of the formation of the solar system) due to different treatments of infall is only about 15%, much less than that caused by different IMF choices.

The different behaviors of nitrogen in our models are illustrated in Figure 7 (a). We focus on N behavior at the same low metallicities which are observed extragalactic H II regions, and we include data from Pagel et al. (1992) and Skillman et al. (1994). While there is a large scatter in the data, the trend seems better fit by the models with mixed and the purely primary N. Note in particular that the mixed model does a surprisingly good job at the low metallicities. While the fit is to somewhat an accident of our particular choices for the primary/secondary N mix, it is amusing that the lowest metallicity region—which dominates the $Y_p$fit—comes out as well as it does.

The effect of secondary N on the $Y-$N relation is shown in Figure 7 (b), with the $Y-$O relation in Figure 7 (c). For these plots we have chosen models with elevated ^{4}He production, as this provides a better fit to the data, (§4.4.). The models with pure secondary N do indeed have a sharp increase at low metallicities; this arises from helium production in concert with suppressed N production due to the dearth of its seed nucleus C. Note, however, that the low metallicity $Y-$N relation is much more gentle for the case of partial primary and partial secondary N, the case that best fits the observations. The small departure from linearity in this case suggests that linear $Y-$N fits might not be in much error (§4.4.).

4.2. Global Constraints and Model Uncertainties

Having sketched some of the effects of the global model features, we now explore the range of abundances calculated for the initial suite of candidate global models and for the subsequent set of light element models. In finding the range of light element predictions we obtain a rough quantitative estimate for the model uncertainties inherent in the chemical evolution arguments used to deduce the primordial abundances of these elements.

We created the suite of candidate global models by making every allowed combination of features shown in table 2; this produces 1184 models. We ran all of these models, using one particular choice of light element model features. Results of this run are summarized in Figure 8 (a), in which we plot the abundances for each element, calculated at the birth of the solar system ($t_\odot$, 4.6 Gyr ago). Abundances are expressed in terms of the logarithm with respect to the observed solar abundance A/H$_\odot^{\rm obs}$:

$$[A/H] = \log \frac{A/H(t_\odot)}{A/H_\odot^{\rm obs}} \qquad (31)$$

(solar data is from Geiss (1993) for D and ^{3}He, and from Anders & Grevesse (1989) for all other elements). Note the wide range of all calculated abundances, including those for the light elements. Indeed, the only element without large excursions is ^{4}He, since it is the only element

whose solar system abundance is mostly primordial and so on this scale does not show large effects of uncertainty due to chemical evolution.

We now impose the constraint that our models reproduce the solar abundances of CNOFe within ± 0.4 dex, i.e. within a factor ~ 2.5. While this may seem generous, recall that we have not tuned our models to fit any elements in particular, yet even models that do so (e.g. Timmes, Woosley, & Weaver 1995) allow themselves a factor of 2. Also, the scatter in the age–metallicity relation suggests that there is a large spread of metallicities at a given time. We must be aware of the possibility that the solar data will not necessarily reflect the average at the solar neighborhood, so we do not want to exclude potentially good models on the basis of a single element.

When we impose this constraint on the candidate global models, we reduce the field to 267 models. A plot of the solar abundances for these models appears in Figure 8 (b). The light element uncertainties are seen to be greatly reduced, while still being considerable. The spread in D and ^{3}He is of particular interest, as we have not added any model variance in the D or ^{3}He evolution (e.g. different η or different low mass ^{3}He stellar processing). Thus the range abundances of D and ^{3}He reflect only the variance due to the global models. Quantitatively, we have a variance in D of a factor of 2.0. As we have noted, this is very close to the range due solely to the IMF, which we see to be the largest source of D variation. We also find ^{3}He to vary by a factor of 1.9, and ΔY to vary from 0.018 to 0.046.

As discussed in §3.1.4., we also constrain the models with infall to fit the G-dwarf distribution, again with a loose tolerance. We focus on the ability the models to avoid a large number of low metallicity G-dwarfs, comparing our results to the data of Rana (1991). A plot of the distributions for one model that passes and one that fails can be found in Figure 9. As discussed in §3.1.4., we do not apply this constraint to models without infall. With the constraint, we reduce

the number of models to 114, with 28 of these being infall models.[5]

We now have a set of 114 chemical models which fit the solar system behavior within the tolerances we have set. To investigate the range of light element evolution allowed by these models, we run each of them for 18 different varieties of light element evolution models. Note that this is far less than the possible number of allowed combinations of options for light element evolution. However, to run such a number of models would not only be tremendously time consuming, but it would also be redundant. We do not expect, for example, the various LiBeB yields in the cosmic ray models to depend on η in an important way compared to the other large uncertainties. Thus we have reduced the number of light element variations to only those 18 most likely to give interesting results. In total, therefore, we have 18 light element models $\times$ 114 global models = 2052 overall models to run.

Results for the 2052 models appear in Figure 10. Notice that a problem has developed in the CNO abundances—some light element evolution features have affected them. The problem arises due to the extra CNO contribution from massive stars with mass loss. For consistency with our protocol of constraints, we again constrain the models to give CNOFe within ± 0.4 dex of solar.

This reconstraint reduces the number of models by about 25% to 1460. Results appear in Figure 11 (a). Note that even for this smaller number of models, the range in the light element abundances remains very similar, although with fewer models the extreme variations are less populated. The models depicted in Figure 11 (a) will

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[5] The only single features clearly important in determining whether a given model would pass or fail were the IMF slope and mass limits. We found no models with $x = 2$ survived; for $x = 2.35$, no models with the widest mass limits passed; for $x = 2.7$, mostly the models with the smallest mass limits passed. Aside from the IMF, no other model feature was prominent in either its presence or absence from the passing models.

be the basis for the analysis in the sections below, except where otherwise noted; we will refer to these as the "final set" of models.

The source of the gap in the D abundances in Figure 11 (a) becomes clear when we plot the results for a particular η. As Figure 11 (b) shows, the D gap arises due to the very high initial D at $\eta = 2 \times 10^{-10}$. It is also clear that it is hard for our models to reduce such a high initial D to its presolar abundance, a point we will explore in more detail in the next section. Note that ^{3}He also shows sensitivity to its initial abundance.

The spread in the light element abundances, particularly the spread beyond that in the "control" group CNOFe, gives a sense of the spread due to model dependence. Note, however, that one must be careful in interpreting these limits. While we have tried a broad sample of models, the lack of a firm theory of star formation leaves open the possibility that one might construct other ones. Moreover, recall that we have not used chemical evolution to get limits on the primordial abundances. Instead, we have assumed BBN yields and seen how well these fit the solar system. What one really would like to know is not this but rather the opposite case. Namely, one would construct models with variable primordial light element abundances (i.e. not necessarily given by BBN values at a given η), and constrain these models to reproduce the solar system values for the light elements and everything else. The resulting spread in the primordial abundances would be an indication of the model uncertainties incurred when setting limits on η. This procedure suffers from an exclusive reliance on the veracity of chemical evolution models. Furthermore, we may estimate the model ranges such a calculation would give, as we now discuss.

Table 3 summarizes the model variations in the light element production for each η. We see that, for all η, D varies almost exactly by a factor of 2 and ΔY varies by 0.063 units. The variation in D+^{3}He depends somewhat on η, but

is small, varying by $< 20\%$ from ~ 2.8. The η-independence of these quantities are evidence that the range model variations does not depend on the light element initial values. This gives us confidence that, were we to do a light element evolution study by allowing the primordial values to float (as described in the preceding paragraph), we would find similar ranges. That is, we expect the reduced quantities in table 3, namely D_{max}/D_{min}, $(D+^3He)_{max}/(D+^3He)_{min}$, and $Y_{max} - Y_{min}$, to be generic estimates for the chemical evolution model ranges of these elements. Of course other selections of model features will give different numbers, but these results at least give a quantitative estimate for the model dependence within a large scale of models, and serve as a benchmark for others.

We emphasize that the D variation is due solely to variation of global model features, in particular the IMF and also infall. Thus to reduce the variation in D requires better chemical evolution modeling as a whole. By contrast, the span in the ^{3}He and ^{4}He ranges are due in part to the effects of different stellar processing features for these elements. Improvements in these features could reduce the ranges given, independently of improvements in the global chemical evolution framework.

It is encouraging to note that the variations in D, ^{3}He, and ^{4}He, while large, are significantly smaller than the 0.8 dex (i.e. factor of 6.3) overall variations allowed for CNOFe in our global model constraints. This implies that the light elements through ^{4}He have less model uncertainty than do the heavy elements. However, the relationship between the heavy and light element uncertainties is not necessarily a simple one.

To get a feel for the connection between the heavy and light element model variations, we examine the effect of tightening the fiducial tolerances in the heavy elements. Specifically, we now require that our models reproduce solar CNOFe to within ± 0.3 dex, i.e. to within a factor of 2. Results appear in Figure 12. The reduction

in the number of models is drastic, going from 1460 models at ± 0.4 dex tolerance to 479 models at ± 0.3 dex tolerance, a 66% reduction in models from only a 20% tightening of the constraint. One can understand this large reduction by noting that our heavy element yields generically are scattered around the solar levels. For example ^{12}C is produced on "average" above the solar level, while ^{16}O is produced in a large range centered below solar; in other words, the O/C ratio is typically subsolar. Thus our models have some range of error in reproducing O/C$_\odot$, and will never have all abundances any more accurate than this range. Indeed, demanding simultaneous fitting of both abundances for progressively smaller tolerances would eventually eliminate all models when the constraints become too stringent given the heavy element yields we have used.

Despite the large reduction in number of models which survive the stronger heavy element constraint, the spreads in the light element abundances remain roughly the same. This reinforces the notion that the CNOFe spreads are to some degree decoupled with the light element spreads. More quantitatively, we find for these models that D varies by about a factor of 1.76 and ^{3}He by a factor of 3.1, values closer to the range for models with the ± 0.4 dex CNOFe tolerance than the 20% tightening would naively allow. In other words, by making the variation in CNOFe smaller by 20%, we do not reduce the variation in D and ^{3}He by as large a factor. Thus the model variation in the heavy and the light elements are not related by a simple linear scaling, but instead the relationship between the two seems to be more subtle.

4.3. Correlations in Light Element Evolution

We now study in detail the results from the set of 1460 models comprising the "final set" having all allowed light element variation while also fitting CNOFe. In particular, we will be interested looking at D, ^{3}He, and ^{4}He (pre)solar and

ISM predictions simultaneously to see better the correlations in their evolution and to show the interplay of the observational constraints.

In Figure 13 we plot the presolar D abundance versus the presolar ^{3}He abundance for each model. One notices immediately that the choice of η (different shaped points) is the largest effect controlling the presolar abundance of both elements. Furthermore, the correlation between D and ^{3}He is evident: the negative slope in the lines traced out in the Figure points out the well-known tradeoff that D destruction leads to ^{3}He production. The steepness of these lines (which radiate out from the primordial values at each η), depends on the degree of low mass destruction or production of ^{3}He. The highest of the three "prongs" for each η comes from the case of ^{3}He production in low mass stars. The middle prong is the "break even" case, and the lowest one from net ^{3}He destruction. We clearly see that the ^{3}He presolar yields are very sensitive to the low mass stellar behavior, which we have only varied by 30% either way from break even. Our models thus confirm the results of Vangioni-Flam, Olive, & Prantzos (1994) and Olive et al. (1994), and extend them to include a very wide range of models.

Figure 13 also includes lines marking the $2-\sigma$ variation in the observed presolar D and ^{3}He abundances, as calculated by Geiss (1993). The Figure provocatively shows that none of the low-η, i.e. high initial D and ^{3}He, models are able to fit the presolar data to within this tolerance. Indeed, the majority of the moderate-η models also seem to fail. At face value, this might suggest that low values of η are ruled out and that higher ones are favored. The question is whether this strong conclusion holds up under scrutiny.

As we have discussed in the previous section, it is a subtle matter to use our models to constrain the possible BBN abundances. We have not included all possible chemical evolution models (e.g. IMF and SFR schemes) and one cannot conceivably do this, given the freedom one has

to adjust each. Also, recall the observed scatter in Fe abundances for stars which are presumably coeval. If the presolar D is a low variation to the mean we have calculated, we would be incorrectly constraining our results. Finally, as we have noted, significant uncertainties and assumptions underlie the entire chemical evolution framework we have adopted (and variants thereof). Even if one could show that if no such models work to explain the low-η D and ^{3}He, such a conclusion would not be founded upon a first principles calculation, since none exists.

Nevertheless, it is not terribly surprising that $\eta = 2 \times 10^{-10}$ fails for our study. Indeed, as we have noted, this value is outside of the "sensible" range in η given by Copi, Schramm, & Turner (1995), although it is within the "extreme" range. Also, while chemical evolution models are uncertain, as we have argued, the D evolution is perhaps the most certain of any nuclide considered. Thus there is no obvious reason to disbelieve the D results, particularly when other abundances can be fit reasonably well.

Furthermore, while chemical evolution is known to be uncertain, our results as presented help one to quantify the size of the model dependence. At $\eta = 2 \times 10^{-10}$, the D abundance varies by a factor of ~ 2; but the lowest D abundance for this η is still a factor of ~ 1.5 away from upper limit to the solar system observation (and with an undesirable ^{3}He abundance at that). It is incorrect to think of the spread in our results as some sort of distribution around a mean model, as we have not just performed smooth parameter variations, but we have also changed whole prescriptions for, e.g., the SFR. Thus it is hard to say more than that the presolar value of D for BBN with $\eta = 2 \times 10^{-10}$ is not reachable in our models and seems hard to reach by models similar to ours.

While presolar D is a strong constraint on models with high η, we see that presolar ^{3}He is a very strong constraint for all η. Because ^{3}He survives processing of both D and ^{3}He, its abun-

dances grows with time.[6] However, the presolar ^{3}He is not much larger than its primordial level, and so demands that models produce very little ^{3}He. We see that models with high η, i.e. with the lowest initial ^{3}He, are the only ones that can fit the presolar constraint. Within these, the different ^{3}He processing becomes important, with ^{3}He production in low mass stars often overproducing ^{3}He. Since this low mass behavior nevertheless seems to be demanded by planetary nebula data, the data seems to favor models which have very little processing in general.

We may also examine correlations between the calculated (pre)solar abundances of other pairs of light elements. Figure 14 shows D and ^{4}He (pre)solar abundances. Here again, the largest segregation of points is due to the different initial D abundances for different η. The shapes of the point distributions are similar for the different η values, though the higher η regions are more compressed. Note that the regions are shifted upwards for increasing η, reflecting the higher $Y_{\rm p}$ for these models. However, it is clear that the shift due to different $Y_{\rm p}$ values is a much smaller effect than the spread due to chemical evolution effects, which give ΔY anywhere from 0.04 to 0.08.

For a given η, the D–^{4}He correlations are not as immediately clear as for the D–^{3}He plot. One can understand the trends by separating the points according to the ^{4}He model features to be varied, namely the $8-12$ $M_\odot$ yields and the presence of mass loss in massive stars. For models with no extra ^{4}He yields (i.e. using just the adopted yields without mass loss and interpolating between $8-10$ $M_\odot$), the predictions appear as the lowest band of points; data for these models alone appears in Figure 14 (a). These models do indeed show an appreciable correlation, as we have argued they should: D depletion is the measure par excellence of the amount of material pro-

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[6]This is true for sufficiently early epochs and for sufficiently high g_3, both of which are the case in our study.

cessed through stars, and ^{4}He production is one consequence of this processing. Thus we expect and find a negative correlation for the "standard" case.

Figure 14 (b) shows the models with one or both of the helium enhancing features. We see here, like the D-^{3}He plot, there are three "prongs" for each η. These correspond (for the most part), to the three different cases of "nonstandard" ^{4}He yields. The lowest trend is for mass loss only, the intermediate (and some of the highest) prong is due to high $8 - 10$ $M_\odot$ yields, and the highest points come from a combination of the two. Here again, we see the negative correlation we expect. We also note that many models appear to fit the solar ^{4}He. Indeed, while some are high and some low, a good number of both standard and nonstandard ^{4}He yields give a successful fit.

We should emphasize that the physics that underlies the D–^{4}He correlation is different from that behind the D–^{3}He correlation. In the latter case, we have demanded that D be made into ^{3}He, and so we have explicitly built in the correlation between the two. The relation between D and ^{4}He is somewhat less direct, as Galactic ^{4}He does not come from primordial D, but rather the destruction of D happens in processes which also produce ^{4}He independently of the level of D. Finally, having found correlations between D and ^{3}He as well as D and ^{4}He, it one also expects correlations between ^{3}He and ^{4}He production. These exist but are more complicated at there is an interplay of both the ^{3}He and ^{4}He evolutionary features. The basic trend, however, is unsurprising: a correlation between the two.

Having examined the correlations between different model calculations at the epoch of solar birth, we now turn to correlations at the present epoch. Here the underlying physics is the same and so we of course expect similar relationships between the model yields, but the observational constraints are different, and in the case of D, more severe. In Figure 15 we plot D versus ^{3}He

for the present epoch, i.e. we plot the abundances calculated for the ISM now. One notices the same features as the plot for the solar system, now smeared out with the intervening evolution. For ^{3}He the observational situation is uncertain as the data shows a large dispersion; if one included the range of abundances spanned by the observations it would not constrain any of these points. This contrasts with the power of the presolar ^{3}He, which offers a tighter constraint than the presolar D.

The D observations in Figure 15 are from the *HST* observation of Linsky et al. (1993), with the $2 - \sigma$ variation shown. (Unfortunately, we are unaware of accurate ^{4}He abundances in the ISM, and so we are unable to constrain our models with this isotope.) If this number is to be taken seriously, it is a strong constraint on chemical evolution models. Here again, we find that the low-η points are very disfavored by the D observations, with the lowest points in this regime needing to move at least a factor of 2 to meet the data. On the other hand, the spread in these points is larger than the presolar data, being almost a factor of 3. Recalling the caveats above, we note that our models have trouble fitting the ISM D abundance given an initial D from a low-η BBN model.

The stringency in the observed presolar and ISM abundances for D, illustrated in figures 13 and 15, bears further investigation. As we have noted, D is the nuclide whose chemical evolution is the simplest, and thus we can expect a given model to calculate D evolution the most accurately of all the nuclides. It is thus interesting to ask what models survive the constraint that both the presolar and ISM D observations are fit to within $2 - \sigma$. As is clear from figures 13 and 15, the more stringent of the two cuts is comes from fitting the ISM observation.

Demanding that models fit presolar and ISM D reduces the number of models from 1460 to 231. In Figure 16 we plot the calculated presolar D versus ^{3}He for these models. Comparing these

results to those of Figure 13, we see that some of the models which could otherwise fit the (also stringent) presolar ^{3}He are removed. However, many of these models still remain. Encouragingly, most of the models which successfully fit the presolar and ISM D also fit the (pre)solar ^{3}He and ^{4}He.

While it is not our purpose to find the "best models" for light element chemical evolution, it is useful to consider the reason for the success of the models which fit the observed (pre)solar D, ^{3}He, and ^{4}He abundances, as well as the ISM D. Most importantly, these models all have a high η and so begin with a very small amount of D and ^{3}He—indeed, at D/H$_p$ = 2.3×10^{-5}, they are already smaller than the $2 - \sigma$ upper limit to the presolar abundance. Thus these demand very little stellar processing of D to reduce it to the observed abundance, and therefore need to avoid producing much ^{3}He.

Interestingly, these models have very few other strong identifying characteristics. All low mass ^{3}He prescription are represented in an even way, as are all combinations of ^{4}He prescriptions. The global characteristics include all SFR prescriptions and infall as well as closed box models. The two steepest IMF slopes are both represented as well, although the only mass limit range is the most restricted one (type 2 in table 2).

We are left to conclude that D evolution is a strong constraint on chemical evolution models, but also that models meeting this constraint can still fit ^{3}He and ^{4}He. We have demonstrated that comprehensive evolution of the light elements is possible within our adopted framework of chemical evolution, although this framework allows for substantial variation in its results according to the adopted model prescriptions and parameters.

4.4. Additional Considerations

4.4.1. D

The program of observation of D in quasar absorption line systems, now just beginning, is likely to revolutionize how BBN is observationally constrained. The power and potential importance and of this method impel a close scrutiny of its assumptions. In particular, it is implicitly assumed that the observed D abundance in high redshift systems is a direct record of the primordial abundance. However, by a redshift of $z \sim 3$, the universe it at least 1 Gyr old, and as much as 3 Gyr old (the range comes about since the $t(z)$ relation depends on Ω_0). Potentially, in this amount of time there might be some depletion of D in the protogalaxies that comprise the absorption line systems; the question is how large such depletion could be. Also, it is noted that typical absorption line systems are not as metal poor as that studied by Songalia et al. (1994), Carswell et al. (1994), and Tytler & Fan (1994); it is thus desirable to understand the evolution of D as a function of metallicity.

Regarding the depletion of D in the absorption line systems, we note first that if these are protogalaxies, we would expect them to take some time to form; however, we model the Galaxy once it has formed. Thus an object at a universal age of 1 Gyr would correspond to a time earlier than 1 Gyr on our models (indeed, Timmes, Lauroesch, & Truran (1995) argue for a significant delay before the onset of galactic nucleosythesis). And furthermore, a glance at Figure 5 suggests that the D depletion at very early times is minimal.

To address this question more systematically, we have considered the D depletion in all of our models for universes at a redshift of $z = 3$, where we have made the very conservative assumption that the protogalaxies form immediately and so time in our models corresponds to universal time. The correspondence between the observable z and the time t in our models depends on the adopted cosmology, in particular on the value of Ω. To be conservative we have chosen $\Omega = 0$ (i.e. a curvature dominated universe, which has the largest age at $z = 3$), with a Hubble constant of $H_0 = 75$ km/s/Mpc. For each of these cases we run our models to the time corresponding to

$z = 3$ and find the D abundance. To get a sense of how this correlates with the metal production, we plot the D abundance for these models against [Fe/H]. Results appear in Figure 17. The Figure confirms that to a high accuracy ($\gtrsim 20\%$ and usually better than 5%) D may be considered undepleted in these systems. Furthermore, Figure 18 illustrates that for metal poor systems with [Fe/H] $\lesssim -1$, and not necessarily at $z = 3$, the D depletion is minimal regardless of the IMF and infall choices.

4.4.2. ^4He

The results from the previous section make clear that our ^{4}He options can produce very different ^{4}He abundances at the birth of the solar system and in the ISM. We now turn to the issue of how these features affect ^{4}He at low metallicities, in particular how the model features affect the $Y-$CNO relation studied in extragalactic H II regions. We show this relation for all four ^{4}He model options (all at $\eta = 4\times10^{-10}$) in Figure 19. We see that the $8-10M_\odot$ enhancement is effective at increasing the ^{4}He slope, making an enhancement of about a factor of two over the standard case. On the other hand, mass loss is unimportant in this regime. This result is not surprising when one considers that mass loss—at least as we have modeled it following Maeder (1992)—not only increases ^{4}He but also CNO. Thus while the ^{4}He production is enhanced, the tracers are as well and so the net effect on $Y-$CNO is small. A note of caution, however, is that the mass loss calculations are difficult and so the results we use are subject to large uncertainty; also, they are only available for two metallicities. Improved modeling is crucial to help address the issue of the $Y-$CNO relation.

Figure 7 (a) shows effect of secondary N evolution on the low metallicity $Y-$N relation; this relation needs to be well understood to properly extrapolate from the observations to derive the primordial ^{4}He abundance. Oftentimes a linear relation is assumed; we now estimate the error

in assuming a linear $Y-$N relation for different cases of N evolution. For each model, we will try to reproduce the empirical fitting procedure by finding the slope of the $Y-$N curve at a point and then extrapolating to get the estimated primordial ^{4}He abundance Y_p^{est}. We can then compare this to the actual Y_p value in the model and thus compute the error $\delta Y_p = Y_p - Y_p^{\text{est}}$ in this extrapolation procedure.

A potential complication to this procedure is that those models lacking features which enhance ^{4}He (mass loss and enhanced $8-12M_\odot$ production) will have smaller $Y-$N slopes and so will be poor approximations to the extragalactic H II region data. Thus we will also keep track of the average slope one would assume for these regions if one did a linear fit; only the results with a sufficiently high slope will be admissible for this comparison.

We perform the extrapolation at N/H = 50×10^{-7}, i.e. at a region surrounded by data and not at low enough metallicity to itself betray the low-N dropoff in ^{4}He. The average slope is not the one used in the extrapolation but is taken at N/H = 100×10^{-7}. This quantity serves only as a diagnostic to indicate the goodness of fit of the given model's $Y-$N relation to the observed slope. As such, a central point better quantifies the average behavior over the whole region.

In Figure 20 we plot the error δY_p in the ^{4}He linear extrapolation versus the average slope dY/dN one would assume for the region. Note the distribution in slopes; we will only consider models with slopes above 10×10^4, which is a rough lower limit to the slopes derived from the H II region observations. For the models with large slopes, one can immediately discern the trends seen in Figure 7, here writ large. Namely, the models with no secondary N are well fit to a line, showing little error in the extrapolation. Indeed, the errors are fairly evenly spread around zero, with a width of about 0.001 units in the mass fraction. The models with some secondary and some primary N give $\delta Y \sim 0.002 - 0.004$,

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while models with purely secondary N can have corrections as large as 0.015.

Thus the models with exclusively secondary N give $\delta Y_{\rm p}$ at the level claimed by Fuller, Boyd, & Kalen (1991), Mathews, Boyd, & Fuller (1993), and Balbes, Boyd, & Mathews (1993). However, as pointed out by Olive, Steigman, & Walker (1991), the N–O relation in extragalactic H II regions demands some degree of primary N; indeed, Figure 7 fits the data well with a mixture of primary and secondary N. For this case, the error in $Y_{\rm p}$ extrapolation is small. Indeed, differences at this level are unimportant compared to the $\delta Y \simeq 0.008$ observational uncertainty.

It is amusing to note that linear extrapolations of H II region data often give slightly larger values when using N as a tracer than when using O, with the difference being around 0.002 to 0.003 units, just the level of the effect we see in the primary + secondary N models. Clearly more work on this issue is needed, and the good agreement of our mixed N model may be somewhat fortuitous, as we have not explicitly built models of the irregular galaxies in which these observations are made. Nevertheless, we our models suggest that the the effect of secondary N on ^{4}He extrapolation may not be as pronounced as originally suggested.

5.   Conclusions

Our results on the sources of our model sensitivity confirm those of Tosi (1988), as we find that the elemental abundances of our models are most sensitive to the shape and mass range of the IMF. Indeed, the models we tried were incompatible with a very shallow IMF slope for any mass range; this was the only model feature that was completely excluded by our solar system constraints. Furthermore, we find that our "best" models, those fitting the solar CNOFe as well as solar and ISM observations of D, select an intermediate IMF slope, namely the Salpeter (1955) value. We also find significant sensitivity to the

presence and timescale of infall. However, infall itself is strongly constrained by the low metallicity G-dwarf distribution. Finally, we do not find our results to depend strongly on the choice of the star formation rate.

We also follow Tosi (1988) in finding that our results are very sensitive to the stellar nucleosynthesis yields employed. We emphasize in particular the effect of systematic effects such as mass loss which are not included in many models. In our models mass loss can be an important source of ^{4}He as well as CNO; more detailed calculations of this affect are needed to address the important question of the ^{4}He contribution of high mass stars. We also find that our results are sensitive to the behavior of stars in the poorly understood $8 - 12\ M_\odot$ range; this too provides an important uncertainty for ^{4}He production. Work on these objects would be as useful as it will be difficult. Finally, we stress the need for a good understanding of the ^{3}He yields in stars of all masses. While it would be quite useful to update the Iben & Truran (1978) calculation of ^{3}He production in low mass stars, it is as important to understand ^{3}He behavior in high mass stars, with emphasis on the effect of winds.

Regarding the viability of light element chemical evolution, it is an important point simply that we do find models which can fit solar and ISM constraints on D, ^{3}He and ^{4}He. This important "existence proof" demonstrates that the framework we have adopted, despite its roughness, is able to fit the data. Thus a consistent picture of light element evolution, from the early universe to the present day, can be drawn.

From the perspective of BBN, it is reassuring that we find the effect of different η so important for D and ^{3}He chemical evolution. We see that the global chemical evolutionary models we have examined, while having a significant degree of model dependence, nevertheless preserve information about the primordial abundances. Thus we are encouraged in our use of solar and ISM data as meaningful constraints on BBN, with the

caveat that there remain significant uncertainties arising from chemical evolution. Indeed, as we have shown, for D and ^{3}He the model dependences are larger than the observational uncertainties (for solar D and ^{3}He as well as ISM D). Thus BBN constraints coming from the solar and ISM D and ^{3}He are useful but should be regarded with caution.

On the other hand, we do not find D to be subject to significant depletion or chemical evolutionary uncertainty at early epochs probed by measurements of quasar absorption line systems. We find the D abundances at this epoch to be a clean indication of the initial D abundance (and thus of η); this result is independent of the particular cosmology chosen. Thus more observations of these systems[7] should eventually be able to put strong constraints on η and thus on $\Omega_{\rm B}$.

In contrast to the case of D and ^{3}He, we find ^{4}He evolution to be dominantly sensitive not to its initial abundance but to the uncertainties in its stellar production. However, we do find that within these uncertainties it is possible to find ^{4}He evolution at low metallicity which can fit the $\Delta Y/\Delta$N and $\Delta Y/\Delta$O slopes better than standard models would indicate. Further, we find that the low metallicity $Y-$N trend is indeed sensitive to the prescence of secondary N. However, our models which best fit H II region N and O data require some degree of primary N and consequently do not lead to large deviations from linear $Y-$N relations. This issue, however, merits further study.

Future work in light element chemical evolution in part would refine the framework adopted here, most importantly by the adoption of different models of the halo phase. One might eventually hope to create a self-consistent stellar and chemical evolution scheme for the light elements, along the lines of the approach of Timmes, Woosley, & Weaver (1995). However, there is ul-

timately a need for a more fundamental understanding of star formation and its relation to the dynamics and chemistry of the Galaxy. In particular, an understanding of the physics behind the SFR and the IMF would be of tremendous value. Nevertheless, while uncertainties in chemical evolution remain, it is encouraging that they still allow models for the light elements to teach us about about BBN.

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[7] Note however that deriving D abudances from the observations may not be simple (Levshakov & Takahara 1995)

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Table 1: Galactic Sources for the Light Elements

NUCLIDE	PRODUCTION SITE	CONTRIBUTION TO OBSERVED ABUNDANCE
D	none	not applicable
^{3}He	low mass ($\lesssim 3M_\odot$) stars (?)	$\gtrsim 50\%$ (?)*
^{4}He	stars of all masses	$\sim 10 - 20\%$

* ^{3}He abundances in the ISM vary with mass of
H II region

Table 2:   Global Model Features

Model Feature	#	Descrption
SFR	1	$\nu\sigma_{\mathrm{tot}}$
	2	$\nu\sigma_{\mathrm{g}}$
	3	$\nu\sigma_{\mathrm{tot}}\mu^2$
	4	$a\exp(-t/\tau)$, $\tau = 7.5$ Gyr
	5	$\tau = 15$ Gyr
IMF shape	1	$\phi \propto m^{-(1+x)}$
	2	from PDMF (Scalo 1986)*
IMF slope	1	$x = 1.0$
	2	$x = 1.35$
	3	$x = 1.7$
IMF mass limits	1	$(m_l, m_u) = (0.2,100)$
	2	$(m_l, m_u) = (0.1,60)$
	3	$(m_l, m_u) = (0.4,30)$
infall		$f = f_0\exp(-t/\tau_{\mathrm{inf}})$
	1	$f_0 = 0$
	2	$\tau_{\mathrm{inf}} = 2$ Gyr; 99.9% σ_{tot}
	3	$\tau_{\mathrm{inf}} = 4$ Gyr; 50% σ_{tot}
	4	$\tau_{\mathrm{inf}} = 4$ Gyr; 99.9% σ_{tot}
N yields		$X^{\mathrm{ej}} = X_0^{\mathrm{ej}}(\alpha + \beta(X_C/X_C^{\odot}))$
	1	$\alpha = 1$, $\beta = 0$
	2	$\alpha = 0.5$, $\beta = 1$
	3	$\alpha = 0$, $\beta = 1$
	4	$\alpha = 0$, $\beta = 2$
SN Ia normalization	1	$\lambda = 0.05$
	2	$\lambda = 0.007$

*Only calculable for exponential SFR

Table 3: Model Ranges for Solar D, ^{3}He, and ^{4}He Production

η	10^5 D_{min}	10^5 D_{max}	10^5 $^3He_{min}$	10^5 $^3He_{max}$	Y_{min}	Y_{max}
2×10^{-10}	5.90	11.9	4.26	18.9	0.250	0.313
4×10^{-10}	1.82	3.66	2.00	7.25	0.258	0.322
6×10^{-10}	0.912	1.83	1.44	4.26	0.262	0.326

η	D_{max}/D_{min}	$(D+^3He)_{max}/(D+^3He)_{min}$	$Y_{max} - Y_{min}$
2×10^{-10}	2.02	3.03	0.063
4×10^{-10}	2.01	2.86	0.064
6×10^{-10}	2.01	2.58	0.064

FIGURE CAPTIONS

1. Big bang nucleosynthesis yields as a function of the baryon-to-photon ratio η. The ^{4}He abudance is plotted as a mass fraction Y_{p}; all other elements given as number relative to hydrogen. Open points show abundance used for the three values of η considered in this study. The dashed lines give the bounds for the Copi, Schramm & Turner (1992) "sensible" bounds on η.

2. (a)The different adopted star formation rates, plotted as a function of time for models without infall. (b) The star formation rate $\psi = \nu\sigma_{\mathrm{tot}}$ plotted for the different models of infall. (c) As in (b), for $\psi = \nu\sigma_{\mathrm{tot}}\mu^2$.

3. The (a) total and (b) gas mass surface densities as a function of time. Plotted for $\psi \propto \sigma_{\mathrm{tot}}$ with an IMF slope of $x = 1.35$, with the four different infall prescriptions.

4. Iron abundance as a function of time for different star formation rates. Data are from Edvardsson et al. (1993).

5. For three different IMFs and SFR 5, evolution of (a) D, (b) ^{3}He and D+^{3}He, and (c) ^{4}He.

6. D and ^{3}He evolution for the infall prescriptions. Note the enhancement of D and dilution of ^{3}He in the infall models.

7. Low metallicities abundances of ^{4}He, N, and O. data is for extragalactic H II regions, as reported by Pagel et al. (1992) and Skillman et al. (1994). (a) N versus O. (b) The ^{4}He mass fraction as a function of O/H at low metallicities. (c) ^{4}He as a function of N/H for different treatments of N evolution.

8. (a) Plot of the yields for the full set of 1184 candidate global models. Calculated for $\eta = 3 \times 10^{-10}$ and for a particular set of light element features. Abundances are calculated for the birth of the solar system and compared to observed solar abundances. Solar abundances are from Geiss (1993) for D and ^{3}He, and from Anders & Grevesse (1989) for all other elements; ± 0.4 dex limits to the solar abundances are indicated by the dashed lines.
(b) Yields for the 267 candidate models having CNOFe within 0.4 dex of solar abundances.

9. The G-dwarf distribution in two candidate global models. The model in (a) passes, the one in (b) does not; see discussion in text. Data is from Rana (1991).

10. Solar abundances for the 2052 allowed global models with all light element model variations. No cut has been made for CNO overabundances due to mass loss in massive stars (see discussion in text).

11. (a) As in figure 10, with the constraint that CNOFe be within 0.4 dex of solar; 1460 models displayed. Note that imposing this constraint makes little change in the light element ranges.
(b) As in figure 11, for models with $\eta = 4 \times 10^{-10}$.

12. The data of figure 11 which reproduces CNOFe within 0.3 dex (short dashed line). For comparison, a level of 0.4 dex is shown in a long dashed line.

13. Plot of D and ^{3}He predicted solar abundances for the "final set" of 1460 models.

14. (a) (Pre)solar D versus ^{4}He for models in having no ^{4}He enhancement features.
(b) Solar D and ^{4}He for models having either or both ^{4}He enhancement features.

15. Calculated abundances of D and ^{3}He at the present epoch. As discussed in the text, it is unclear how to use the ISM ^{3}He data as

a constraint; we have omitted it here. We show the D abundance in the ISM due to Linsky et al. (1993).

16. Calculated D and ^{3}He abundances at the solar birth. As in figure 13, but models have been constrained to fit both the presolar and the ISM D abundances.

17. D versus [Fe/H] for our models, plotted at a galactic age equal to universal time at $z = 3$. The redshift-time relation is derived for a cosmology with $\Omega = 0$ and $H_0 = 75$ km/s/Mpc. Dotted lines indicate primordial levels for the three adopted η values.

18. D versus [Fe/H] for the models of figure 5 having different IMFs and for the high infall model 4 of table 2. Regardless of these model features, the D evolution at low metallicity is minimal.

19. Mass fraction of ^{4}He plotted as a function of (a) C, (b) N, and (c) O abundance in the low metallicity regime. Solid line: no ^{4}He enhancement; dashed line: $8-12\ M_\odot$ ^{4}He enhancement; dot-dashed line: mass loss; dotted line, both enhancements. For $\eta = 4 \times 10^{-10}$.

20. The error $\delta Y_{\rm p}$ made in extrapolating the $Y-$N relation linearly to determine the primordial ^{4}He abundance. Plotted for each model as a function of the average slope dY/dN of the $Y-$N relation at low metallicity. Open circles are for models with only primary N, triangles are models with mixed primary and secondary N, squares are models with low secondary N and hexagons are for models with high secondary N. Horizontal line indicates level of observational error; vertical line indicates approximate level of minimum slope for H II region data.